THEORETICAL STUDIES OF THE RS CANUM VENATICORUM STARS

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I. INTRODUCTION AND SUMMARY

The aim of this work has been to improve our theoretical understanding of activity in RS CVn stars. Specifically, we set out to develop models for chromospheric structure, and to examine the role of magnetic fields both in the photosphere as well as in the chromosphere and upper atmosphere. As an extension of the work on RS CVn stars, we also wished to examine the case of the T Tau stars from the same points of view.

In the course of the work done under this grant, we have made significant progress in using the properties of magnetic field loops to unify our understanding of atmospheric structure in RS CVn stars. However, the concepts developed in the case of these stars now appear to be applicable over a much broader region of the HR diagram. The key to our new understanding is the <u>absence</u> of <u>stable</u> magnetic loops in the atmospheres of late-type giant stars. This is a feature for which our experience with the solar atmosphere did <u>not</u> prepare us well. If our new understanding is correct (and we will bring together a broad diversity of observational facts to support it, see below), then the <u>atmospheres of RS CVn active components are qualitatively distinct from the solar atmosphere</u>. This conclusion has an important bearing on research into the broad area of "solar-stellar connections": although detailed knowledge of solar physics and phenomena undoubtedly helps in understanding activity in certain types of stars (e.g. main sequence G,K, and M stars), it will be important to resist the temptation to apply all "solar concepts" universally across the HR diagram.

II. CHROMOSPHERIC MODELS

We have developed a chromospheric modelling code and applied it to calculate flux profiles of H_α , H_β, and H_γ in stars which lie in the region of the HR diagram occupied by RS CVn secondaries. The aim is to undertake first of all a systematic study of Balmer line profiles in order to determine their usefulness as chromospheric diagnostics in these stars. Our work has consisted of three parts: (a) code development for use at Bartol; (b) specific application to λ Andromedae; and (c) computing a grid of chromospheric models which can serve not only to interpret RS CVn Balmer Line profiles, but also in a broader context of understanding the empirical relationship which is known to exist between the width of H_α and stellar luminosity.

(a) Code development

During my stay at Sacramento Peak Observatory in 1978, I had collaborated with Dr. L. E. Cram in developing a non-LTE chromospheric code for a six-level hydrogen atom. At that time, the main interest was to calculate Balmer line profiles in red dwarfs. The code begins by specifying a photospheric model from the literature. The highest point in the published model is taken to be the temperature minimum in the atmosphere, and we superpose on top of that photosphere a rise in temperature upwards into the chromosphere. The temperature is supposed to rise to a value of T_c (typical of the "top" of the chromosphere) at a mass loading m_c. Usually, T_c is taken to be 8000 K, since at that temperature, hydrogen is beginning to ionize appreciably in the atmosphere, and it becomes difficult for the chromosphere

to keep itself cool in the presence of mechanical energy deposition (whose source we do not specify). Above T_c, the temperature is assumed to rise very rapidly, typically reaching 2×10^5K in an interval of only 0.25 in log m. The temperature structure is taken to be linear in log m between m_c and the temperature minimum.

To start the code, a value is chosen for m_c and T_c, with a given photospheric model. An approximate evaluation of the non-LTE departure coefficients for the two lowest levels of the hydrogen atom is then used in order to calculate a preliminary solution of the hydrostatic equilibrium chromosphere. This serves as input for the main program which solves the radiative transfer (R.T.) problem for a hydrogen atom with five bound levels plus continuum. In this code, we solve explicitly for the first three Balmer lines, and for the first three continua of the hydrogen atom. Radiation temperatures are assigned to other transitions and continua, except that the Lyman lines are taken to be in detailed balance. No microturbulence is included: these models are strictly static.

Our solutions for red dwarfs had shown that in such cool stars, essentially no contribution to Hα came from the photosphere. However, this is no longer the case in the stars of interest to us here. In our present work, there are roughly comparable contributions from photosphere and chromosphere, at least in the hotter stars. This requires us to choose our temperature grid in the chromosphere with some care in order to avoid convergence problems. As with the red dwarfs, our code contains an inner iteration (for the RT solution) and an

outer iteration (for hydrostatic equilibrium), because we are dealing with the principal atmospheric constituent (hydrogen), and therefore changes in level population which arise from one converged RT solution to the next can have a serious impact on the hydrostatic equilibrium. (Such an outer iteration does not enter into RT calculations of minor species, such as Mg or Ca.)

Once the main RT convergence has satisfied also hydrostatic equilibrium, we then use the converged model as input to a third program for calculation of detailed flux profiles of the Balmer lines for purposes of comparison with observations. Examples of line profiles from some of our converged models are shown in the Figure 1(a)-(f).

The programs were developed first at Sacramento Peak, and it seemed at first advisable to attempt to convert them to the Delaware machine for use at Bartol. However, all attempts to convert the large RT code failed. I therefore submitted a proposal to Sacramento Peak Observatory requesting time on their PE-3242 machine, which became accessible for remote job entry via a dial-up in June 1981. I was granted about 100 hours of CPU time, and most of the results to be reported here were obtained in this mode of operation.

(b) λ Andromedae: Large Macroturbulence

B. W. Bopp supplied us with four high precision profiles of $H\alpha$ in λ And, taken at various epochs. The observed line profiles are broad, and slightly (few %) asymmetric. This star is especially suitable for study because it is viewed almost pole on to its orbit, and so rotational broadening can safely be excluded when we analyze line profiles.

With T_c=8000 K, we computed several models with log m_c ranging from -∞ to -4. as m_c increases, the absorption feature at Hα deepens and becomes slightly broader at first. But then emission begins to fill in the central regions of the line, and eventually the absorption disappears. Thus, there is a maximum possible broadening of the line in absorption. (We note that it is because Hα responds in this significant manner to changes in m_c that the Hα profiles can serve as a diagnostic of chromospheric conditions in these stars.)

Significantly, all four of Bopp's profiles were much broader (by at least 50%) than we could achieve at maximum absorption strength. Since rotation can be excluded, we tried to broaden the line by macroturbulence. We then found that within ± 2 Å of line center, we could fit the observed line profiles with surprisingly high precision (± 2%), if we used a macroturbulent broadening parameter, ζ, in the range 30-40 km/sec (see Figure 2). (We assumed a Gaussian distribution of macroturbulent velocities.) These velocities seem rather high, and yet the goodness of fit to the observed profiles suggest that they are meaningful.

A different approach to broadening the line is to consider higher pressure chromospheres (i.e. move m_c to deeper layers). However, in this case, in order to prevent Hα from going strongly into emission, we need to reduce T_c to 6500 K. This was the value used by Baliunas et al. (1979) in deriving a previous high pressure model of λ And, and they claimed they could fit the line observed by Kraft et al. (1964) for this star within 10% as regards central intensity and half width. (However, they did not show an illustration of how good the fit

was to the profile as a whole.) There appears to be no physical reason for having T_c=6500 K, unless the mechanical energy deposition rate suddenly increases rapidly above those temperatures. (As mentioned above, 8000 K is more reasonable physically because, with a uniform deposition of mechanical energy, hydrogen will become more than 50% ionized at typically 8000 K.) Nevertheless, for completeness, we also tried a model with T_c=6500 K, and found that indeed, a high pressure model could be calculated which fitted the observed profiles fairly well. (Not as well, however, as before: now the precision of fit was only $\pm$10% (see Figure 2)). Thus, in these models, there is no need to introduce macroturbulent broadening. Our best fits for the λ And profiles turned out to be models with pressures at the top of the chromosphere, p_c, about 0.056 dyn/cm^2 in the low pressure case (including ζ=30-40 km/sec), and p_c=0.4 dyn/cm^2 in the high pressure case.

Can we decide on some other basis between these models? It turned out, after our calculations were finished, that the answer may be yes. Certain spectral lines in the UV have the property that their emission strength ratios are sensitive to densities in the upper chromosphere and transition region (TR). It is true that these line ratios are quite difficult to interpret in terms of absolute densities in the chromospheres, but a differential comparison of the line intensity ratios in two stars ought to be quite reliable as an indicator of density in one of the stars relative to the other. Just such a differential comparison has recently been completed by Ayres et al. (1982) for the stars λ And and α Cen B. For the latter object, a rather reliable model chromosphere already exists, based on intensities

and profiles of many lines in UV and visible portions of the spectrum. From the differential comparison, Ayres et al. concluded that p_c in λ And probably lies in the range 0.05-0.09 dyn/cm^2. Our low pressure model, with large ζ, lies precisely in this range, whereas the high pressure model appears to be excluded (see Mullan and Cram, 1982).

Two other recent studies have appeared in which large macroturbulent velocities (45-100 km/sec) were reported in T Tau stars (Herbig and Soderblom, 1981; Ulrich and Wood, 1981). The activity in RS CVn itself was first described as "T Tau-like" (Hall, 1972) and in fact, there may be a close relationship between T Tau stars and RS CVn stars (see section IV below). Hence, our conclusion of large macroturbulence in the atmosphere of an RS CVn star may be less surprising than we first thought. Moreover, the fact that macroscopically large elements move rapidly through RS CVn atmospheres had been discussed previously by Weiler (1978) in terms of discrete structures ("prominences") in the Mg II k line profiles. The "prominences" in Weiler's data must have been large indeed, since their emission strengths were at times comparable to the integrated emission from the rest of the visible disk of the star. Moreover, the "prominences" which he observed were moving at even higher speeds (up to 250 km/sec) than the speeds which we are proposing here.

So we need to confront the basic question: what source exists for this striking macroturbulence in RS CVn atmospheres? The work which we have been doing on mass loss (see below) suggests one possible explanation. We shall argue that <u>the macroturbulent "elements" are unstable magnetic flux loops</u>. We regard the unifying role played by such unstable flux loops

in interpreting a variety of observational data in late-type giants as the most important conclusion to emerge from the work of this grant.

(c) Grid of chromospheric models: Hα Width-Luminosity Relation (WLR).

For our grid, we used photospheric models as follows: $T(10^3 K)$, log g = 4,4; 4,3; 4,2; 5,3; 5,2; 6,4; 6,3; 6,2. For each photospheric model, we have calculated 4 chromospheric models with T_c and log m_c values as follows: 6500 K, -6; 6500 K, -4; 7500 K, -4; 7500 K, -6. This represents an attempt to span both "high pressure" and "low pressure" models in our grid. Examples of Hα profiles at some of the points in our grid are shown in the Figure 1.

It will now become possible to apply these profiles to interpret observed Hα profiles in RS CVn stars. However, a more immediate question can also be addressed with our results, and this has an important bearing on the overall results of our work, because it stresses once again the importance of velocity fields in late-type stars, and the use of Hα as a diagnostic of chromospheric velocity fields.

The observational relationship between width of Hα and the stellar luminosity was first pointed out by Kraft et al. (1964), and verified by LoPresto (1971). The relationship can be written approximately in the form (where H = half width in Å).

$$\log H \approx -0.047 M_v + \text{const} \tag{1}$$

over a range of M_v brighter than about $M_v \approx 0$ or -2. It is

important to note that the relationship (1) does not extend to less luminous stars, nor to M giants and supergiants, but seems to work only for luminous G and K giants and supergiants. For G and K stars fainter than the above limit, a separate width-luminosity relationship (WLR) seems to hold (Kraft et al., 1964). Part of the difference is that damping wings are beginning to become noticeable in G dwarfs, and this prevents a reliable measurement of the true half-width of the absorption line.

Now, for our models with T_e = 6000 K, log g = 4 corresponds to $M_V \approx +4$ (using Iben's tracks, 1967); while log g = 2 model corresponds to $M_V \approx -3$. Hence, between log g = 4 and 2, equation (1) suggests that H should increase by a factor of about 2.1. Results from our models are shown in the Table.

TABLE 1

H = Half-width of Hα in chromospheric models (T_e=6000 K)

Model	H(log g=4)	H(log=2)	H(2)/H(4)
6500,-6	0.76 Å	0.60 Å	0.8
6500,-4	1.04 Å	0.74 Å	0.7
7500,-4	0.78 Å	0.83 Å	1.1
7500,-6	0.72 Å	0.75 Å	1.0

Our models certainly do not behave in the way the observations indicate: in no case do we find Hα increasing in width by a factor of 2.1 between log g = 4 and log g = 2. It is true that there may be systematic variations of m_c with gravity (Kelch et al., 1978), but these are not sufficient to alter the ratios of H(2)/H(4) to a value as large as 2.1. A similar conclusion emerges from our models with T_e = 4000 K. We cannot explain

the empirical WLR by our static chromospheres, by, e.g. varying the depth of the chromosphere (i.e. m_c). This conclusion is in striking contrast to the explanation which Ayres (1979) has proposed for the WLR in Ca K emission: there, the WLR is due to systematic effects in m_c, rather than in any velocity-related parameter.

Our conclusion therefore seems to be that depth of the chromosphere is not the controlling influence in the Hα WLR. What else can be responsible then, other than velocity fields in the chromosphere? In this regard, the success which we had in using macroturbulence in RS CVn stars comes to mind, and we wonder about applicability of macroturbulence over more general areas of the HR diagram (such as is covered by the sample of Kraft et al.). In this connection, it is worth noting a further important difference between the WLR in Ca K and the WLR in Hα: the former applies essentially in a single-valued way over the entire range of M_v, whereas the latter does not. Kraft et al. point out that the linear relationship in (1) above applies only to brighter stars, and a different relationship apparently holds for fainter stars. The break between the two separate relationships occurs close to $M_v \approx 0$ or -2. Interestingly, values of macroturbulence also show a distinct break at almost the same value of the luminosity (Smith and Dominy, 1979). At lower luminosities, ζ is small for photospheric lines, while for higher luminosities, there is an almost discontinuous jump to larger values of ζ. Now, there seems to be a rather close link between velocity fields in the photospheres and chromospheres of late-type stars (Imhoff, 1977), and so we are led to suspect

that the jump seen by Smith and Dominy in the photospheric velocity fields may be related to the discontinuity in the WLR reported many years previously by Kraft et al., i.e. a discontinuity in a chromospheric velocity parameter.

We are led to the conclusion therefore that Hα profiles in late-type giants may serve as useful diagnostics of chromospheric macro-velocity fields.

III. MAGNETIC ACTIVITY AND MASS LOSS

When we started work on this grant, it had just been discovered from IUE and Einstein data that rapid mass loss in cool giants sets in across a certain boundary line in the HR diagram, which also serves to separate stars with detectable steady X-rays from those with no detectable steady X-rays. During the course of our work in the last two years, it has emerged that this boundary also serves to segregate stars according to other criteria as well. Among these, we wish to draw particular attention to C-13 richness in the atmosphere (indicative of dredge-up processes from deep inside the star), and the behavior of the He 10830 Å line. Thus, the mass-loss boundary which we (Stencel and Mullan, 1980) discovered in the HR diagram, has turned out to be an intensely interesting area for research in several areas of stellar astrophysics.

The feature which had drawn my attention originally to the RS CVn stars was that the mass loss boundary passed through the sub-giants at spectral types K0-K1. Now, it had been known for many years (without any explanation for the fact) that active secondary stars in many RS CVn systems were subgiants of just these spectral classes (Popper, 1970). This suggested

to us a possible connection between rapid mass loss in cool giants and activity. As we shall describe here, this suggestion has turned out to be a fruitful area for research.

(i) Non-thermally driven mass loss

It is important that mass loss becomes rapid among just those giant stars where the X-ray flux (time-averaged) becomes too weak to be detected by Einstein. This forces us to conclude that thermal pressure is not responsible for driving mass loss in late-type stars. Parker's concept of a steady, spherically symmetric wind, with thermal conduction dominating in the energy transport, would lead one to predict that mass-loss rate is very sensitive to coronal temperature: if the temperature drops noticeably (as it appears to do, when one observes giants of later and later spectral types), then, if anything, the mass loss rate ought to be dramatically reduced in going to later spectral types (other things being equal). Precisely the opposite behavior, however, is indicated by the observations. A non-thermal source of the wind must be identified.

During the first month of the grant period, I decided to examine the scenario which is shown in Figure 3. Consider a magnetic loop emerging at a stellar surface (due, e.g. to dynamo activity in the sub-surface gas, which is convective in all late-type giants, and therefore highly conducive to dynamo activity). The way in which such a loop will evolve depends on the properties of both the loop itself (its length, magnetic flux, etc) as well as the properties of the ambient atmosphere into which it emerges. Under certain conditions, it will certainly be possible for such a loop to find a static equilibrium in the atmosphere,

in which case, we are no longer interested in the loop. However, it is also possible for the loop to become unstable, e.g. to a ballooning instability (Low, 1981). In that case, the loop will become greatly distended upwards. Pinching near the base will help to drive magnetic reconnection there, and ultimately a bubble of material will be severed from the star, and ejected as a mass loss "event".

When I presented this concept at the first Cambridge Cool Star workshop, Dr. I. A. Ahmad pointed out that there was already some solar evidence for this type of mass loss process. He and I therefore began a collaborative study of the non-thermal sources of the solar wind.

(ii) Solar wind: magnetic driving?

Skylab data have demonstrated conclusively that the Parker concept for the solar wind is definitely inadequate: the wind is not steady, not spherically symmetric, and certainly is not fastest where the temperatures were highest (coronal holes are cool). Neither can the energy fluxes in the high speed streams be provided by conduction: the conductive energy flux in coronal holes is too small to provide the kinetic energy flux in the streams emerging from the holes. Additional, non-thermal energy is required to drive the solar wind (cf. Brueckner et al., 1977).

Ahmad and Webb (1978) studied X-ray images from Skylab, and examined X-ray bright points (XBP) at the bases of plumes in the solar polar coronal holes. They found that the plumes are the sites of upward flowing material, with substantial mass flux. In fact, taking into account the total number of XBP in the polar

holes (there are typically 5-10 XBP (plumes) per polar hole), they showed that the mass flux was sufficient to explain the entire mass flux in the solar wind. According to Ahmad and Webb, then, the solar wind comes from a small number of discrete sources, any one of which lasts only a relatively short time (less than 1 day, typically). Hence, the solar wind is intrinsically unsteady on a time scale of order one day, and is expected to fluctuate in mass flux by at least 20-30% (assuming a total number of 10-20 "sources" which evolve independently). Statistical confirmation of their hypothesis has been provided recently (during the period of this grant) by Davis (1980), who found a good positive correlation between numbers of X-ray bright points in coronal holes and the mass flux in the solar wind.

This led me to consider the scenario in Figure 3 in more detail in collaboration with Ahmad. We were led to consider a magnetic driver for the mass loss events because (i) XBP are known to be well-correlated with bipolar emerging magnetic flux loops, and (ii) the fact that they are bright in X-rays indicates that energetic processes are already at work in those magnetic loops, presumably related to releasing some magnetic energy in the form of heat. It seemed to us natural to consider tapping some of this released magnetic energy in the form of mass loss. Since reconnection had already been discussed as a possible source of the energy release in XBP (Parker, 1975; Habbal and Withbroe, 1981), and it is well-known that magnetic reconnection leads to ejection of high speed flows (Vasyliunas, 1975), it seemed natural to consider the reconnection process in the evolution shown in Figure 3 as a source of energy for mass ejection. We were able to show that conditions in a loop

are capable of leading to a ballooning instability to start with, and so the upward distending process is consistent with what can happen in the solar atmosphere. We have derived time scales and mass ejection rates, as well as altitudes for dissolution of magnetic bubbles, which were all consistent with solar wind data (Mullan and Ahmad, 1982). At present, the most important aspect of the scenario which requires further research, is just how a loop evolves in a stellar corona subject to a ballooning instability. This is a complicated dynamical problem, requiring numerical MHD modelling. However, it is of sufficiently general interest for the mass loss process from all late-type stars (see below), that it requires detailed attention. The post-doctoral fellow who was hired under this grant (Dr. Wayne Waldron), has been using his plasma physics and astrophysics backgrounds to construct a 2-D MHD code to examine this question.

The major conclusion which has emerged from the work with Ahmad has been the radical re-appraisal to which we have been led for the source(s) of the solar wind. According to this scenario, the sun loses mass in a series of discrete events, each one of which is accompanied by release of sufficient energy to cause local X-ray brightening. Since release of magnetic energy is traditionally considered as the most likely candidate for "magnetic activity" of all kinds, we have here <u>an intimate connection between activity and mass loss</u>. It was precisely such a connection which led us to embark on RS CVn studies, and so the implications of applying what we have learned in the solar case to the stellar case are obvious.

(iii) Solar and stellar coronae: essential distinctions

The key to the scenario for solar mass loss described above is the ballooning instability. It is crucial to note that in the sun, not all magnetic loops are subject to such an instability. Quite the contrary. When a new loop emerges into the solar atmosphere from beneath the surface, it is more likely to enter a long-lived phase of relative stability: closed, quiescent loops are observed to emit X-rays at an essentially unchanging level for days on end in the sun. This is a very important point which our long familiarity with the sun may cause us to overlook. Evidently, the footpoints of a loop are pushed around by the constantly changing convective turbulence near the photosphere, and so in strictness, a magnetic loop evolves according to the MHD equations in time. And yet, most loops can apparently find an almost force-free state, or series of states, through which they evolve in a quasi-steady manner (Low, 1982). At any given time, the loops give the appearance of being stable, because the atmospheric structure is such that they can remain static. It is in this sense that one is led to the basic concept of "building blocks" of the solar corona (Vaiana and Rosner, 1978), i.e. individual loops which remain in existence for long periods of time, and contain essentially all of the coronal emission in X-rays. The term "building block" itself carries a basic connotation of a stable structure. However, when we discuss other stars, it will not be appropriate to carry over the concept of "building blocks" for the magnetic loops, unless these stellar loops can be shown to be stable. The latter is a crucial proviso, as we shall argue below.

In the sun, the fact that coronal loops can be long-lived

and essentially stable, leads to a conclusion which, although obvious, is worth making explicitly in view of what we shall discuss below for other stars: enhanced coronal emission is well correlated with enhanced chromospheric emission. In contrast, if coronal loops are not stable, then there is no a priori reason why chromospheric emission should be positively correlated with coronal emission. In this regard, the following point must be stressed: the boundary in the HR diagram where coronal properties undergo striking alterations (e.g. rapid mass loss, disappearance of steady X-rays), is not accompanied by any sensible alteration in the chromospheric emission strengths of the stars (Mullan, 1981; Mullan and Stencel, 1982). As mentioned above, it is important to prevent solar prejudices from making us overlook the possibility that chromospheric and coronal emission in late-type stars need not be correlated. In fact, just such a lack of correlation has emerged strikingly from He 10830 work by Zirin (1982). The strength of He 10830 absorption is a measure of coronal X-ray flux (because the lower level of the transition is populated via X-ray processes; cf. Zirin, 1975). This strength has been found to be well correlated with the strength of Ca K emission (i.e. chromospheric emission) in G-type stars: our experience with the sun leads us to find this as no surprise. The surprise comes with the K and M stars: for these stars, Zirin finds that the correlation between coronal and chromospheric emissions becomes very poor, essentially non-existent: suggesting that (in Zirin's words) "the late stars of relatively high K-line intensity may indeed have active regions [on their surfaces], but [they] cannot keep the corona from flying away in the stellar wind".

(iv) Helmet streamer stability: solar corona

Pneuman's (1968) work on helmet streamers provides a convenient starting point for analyzing the stability of closed magnetic loops in the solar corona. Helmet streamers are regions of closed field lines near their base, with open field lines surrounding the closed fields. The closed field lines reach out to a maximum radial extent of r_h, while solar wind can flow out along the open field lines, passing through a sonic point at a radial distance of r_s. Pneuman analyzed the steady state case, in which magnetohydrostatic equilibrium obtains in the close field region, and a thermally driven wind exists on the open field lines. Making several simplifying assumptions, in order to keep the discussion analytic, Pneuman derived a simple relation between r_h and r_s:

$$r_h = r_s/2 \qquad (2)$$

I have re-done Pneuman's calculations, relaxing the various simplifying assumptions which he made, in order to determine how (2) is altered from the simplest case. I have found that the relationship in (2) holds over a considerably broader range of conditions than Pneuman assumed, although it is true that the numerical value of the denominator may need to be increased somewhat. Nevertheless, to much better than an order of magnitude, (2) can serve as a useful approximation.

Pneuman found that his steady state situation could exist only as long as the coronal temperature was less than T_{max}. The numerical value of T_{max} turned out to vary (slowly) with an area parameter, β (= ratio of area of closed field lines to area of open field lines). Pneuman had no means of deciding on the

correct value of β, because of his simplifications. Over a broad range in β, (factor of 10^4), he found T_{max} in the range from 0.8 to 3.6×10^6 K. For $T > T_{max}$, (once β is prescribed), no static solution of the helmet streamer exists: all field lines presumably become open. I have tried to make this conclusion more definitive by collaborating with Dr. R. S. Steinolfson (UC, Irvine), who has a 2-D MHD code for the study of coronal transients. In this application, one starts the code with a dipole magnetic field at the center of the sun, and then at t=0, impose a Parker-type solar wind flow. The flow interacts with the field, and vice versa, and eventually approaches a steady state with some field lines closed, some field lines open, and the ratio β is determined by the steady state solution, rather than being a free parameter. For low values of T (coronal temperature), some closed field lines do indeed survive, and equation (2) is satisfied to better than a factor of 2. Moreover, as T increases, a value is eventually reached (between 4 and 4.5×10^6 K), such that at higher temperatures, the steady state solution contains no closed field lines above the solar surface. This is valuable confirmation of the general validity of Pneuman's discussion.

(v) Helmet streamer stability: stellar coronae

In the sun, the only way to force a helmet streamer towards instability is essentially to allow T to increase. However, in other stars, we can also use the gravity to approach instability. That is what I have done in an application of Pneuman's results to the stellar case in general. I used empirical chromospheric pressures as pressures at the base of the stellar coronae (Kelch et al., 1978), and then I assumed the minimum flux coronal concept

(Hearn, 1975). With these, I could ask the question: where in the HR diagram is the condition $r_s = 2r_\star$ satisfied? The results are shown in Figure 4, where I have used Paczynski's (1970) evolutionary tracks to convert masses and radii to points on the HR diagram. The striking result is that $r_s = 2\,r_\star$ along the Mg mass loss boundary, such that stars which are losing mass rapidly have $r_s < 2\,r_\star$, while stars which do not show rapid mass loss, have $r_s > 2\,r_\star$.

The significance of this emerges from considering equation (2). Stars with $r_s > 2r_\star$ have $r_h > r_\star$, just as the sun does, and so closed loops (and steady helmet streamers) may exist in stable conditions in the atmospheres of such stars. Hence, they can be strong in X-rays, low in mass loss rate, and show good correlation between coronal and chromospheric emissions. However, if a star is above the mass loss boundary, it has $r_h < r_\star$, i.e. <u>the last closed field line is inside the star</u>. If dynamo action causes a new magnetic loop to emerge at the surface of the star, then the radial extent of that loop already violates the condition for the last closed field line of a static helmet streamer. Therefore, <u>in such a star, an emerging loop cannot exist in a static helmet streamer type of configuration</u>. Hence, no closed steady loops exist in the atmospheres of such stars. This seems to explain naturally why X-rays from such stars are at a very low level. Furthermore, if closed loops cannot be stable, what can they do when they emerge into the atmosphere? They must evolve dynamically, and this immediately reminds us of the scenario discussed above for mass loss in coronal XBP: we expect that dynamical evolution will enter a ballooning phase, and field loops in late-type giants will be distended upwards in the atmosphere (cf. Fig.3(b),(c)).

In this regard, it is important to remark that greatly distended loops play a major role in interpreting X-ray and IUE data of RS CVn coronae (Walter et al., 1980; Swank et al., 1981: Simon et al., 1980), for entirely independent reasons from the ones which we are discussing here. And as a support of our finding that the onset of extended coronal loops occurs among subgiants at early K type, we may cite recent evidence of X-ray eclipses in the RS CVn system AR Lac, consisting of an early G subgiant and an early K subgiant. The details of the X-ray light curve in the immediate vicinity of both primary and secondary eclipse in this system are such that the X-ray emission around the G star must be confined to a thin corona close to the stellar surface, whereas the K subgiant must have an extended corona, reaching out to radial distances of order 2R(star) (Walter et al., 1981). This system is therefore particularly interesting in the present context since it happens to contain one star on one side of MTTL,(see below) and a second star on the other side of the MTTL, both of which are in a position to eclipse (at least partially) the other.

What eventually happens to a loop which is balloon unstable? So far, a definitive answer cannot be given. However, according to the scenario in Figure 3, magnetic reconnection may ultimately sever the connection with the stellar surface, and a bubble of material will be ejected from the star. In this scenario, then, mass loss among late-type giants is an intrinsically episodic process, with discrete mass ejection events each time a new flux loop emerges at the stellar surface. In this view, <u>the Mg mass loss boundary corresponds to a transition in the overall magnetic topology in the stellar atmosphere from mainly closed to mainly open</u>. Because of that, we call this the "Magnetic Topology Transition Locus" (MTTL).

Above MTTL, mass loss is considered to be driven by magnetic reconnection processes. Since these are generally responsible for magnetic activity also, we propose that magnetic reconnection in unstable loops provide the causal link between activity and mass loss which first drew our attention to RS CVn stars. This is an important point, because if mass loss is being driven by an energetic process of the kind we propose, then <u>transient</u> (non-thermal) X-rays should be observed from time to time in late-type giants. The X-rays may be hard, depending on how strong the reconnecting magnetic fields are (e.g. in RS CVn stars). In fact, rather energetic processes must be occurring in RS CVn atmospheres in order to explain the non-thermal radio outbursts, and the hard X-rays which seem to set the RS CVn systems apart as an especially powerful class of active stars: the hard X-rays from RS CVn systems cannot be easily understood in terms of solar analogies (Garcia et al., 1980). Thus, we make a distinction between quasi-steady (thermal) X-rays which may characterize long-lived loops in solar-like atmospheres, and the intrinsically variable (perhaps hard, probably non-thermal) X-rays which characterize the mass-ejection events in cool giants. In this regard, it is worth noting that the hard X-rays in RS CVn systems are indeed highly variable in time (Swank et al., 1981).

The fact that a source of transient X-rays exists (in our scenario) in cool giants from time to time has an important bearing on the excitation of the He 10830 Å line. This line becomes stronger in absorption, the larger the coronal X-ray flux becomes (Zirin, 1975). In general, therefore, it is expected that in crossing the boundary where mass loss becomes rapid, and (steady) X-rays die away, the He 10830 Å absorption should also die away. In a broad sense, that turns out to be true (Zirin, 1982): He 10830 weakens monotonically in going from G to K to M stars of all luminosity classes from I to IV (dwarfs do not show this behavior), at least when one averages over

broad spectral classes. However, Zirin (1982) has noted that in finer detail, the decline is not monotonic, but shows a peak at K3 among the giants. He o fers no explanation for the fact, but he also points out something which has a bearing in the present context: variability in the profiles (i.e. what Zirin calls "activity") also is most prominent at K3 among the giants. The puzzle of He 10830 excitation in cool giants has been discussed in some detail by Simon et al. (1982), and they also cannot provide an explanation for why the line should be so strong in stars where, on average, the X-ray coronal flux ought to be negligible.

Our scenario provides a natural interpretation: transient X-rays from mass ejection phenomena excite variable He 10830 absorption, and the more pronounced the activity, the more pronounced the He absorption can become. The low photospheric temperatures of late-type giants have the effect that He 10830 absorption is very sensitive to coronal X-ray flux (more so than solar).

(vi) Discrete mass ejection

The He 10830 data also provide evidence which has a bearing on episodic mass ejection. Zirin (1982) has shown that the 10830 line cannot go into emission when seen against a stellar disk without at the same time violating other observational constraints. Therefore, when He 10830 appears in emission (which it does from time to time in some of Zirin's 455 program stars), it must be created in a detached shell of material around the star. Presumably, the shell indicates an episode of mass ejection. As many as 4 discrete mass ejection events can be traced in the records of He 10830 in certain stars (Zirin, 1982), in the course of 10-15 years. Hence, discrete mass ejections seem to be rather common events in stars which have evolved across the MTTL., as our scenario requires.

(vii) Macroturbulence

Unstable magnetic flux loops provide a natural explanation for macroturbulent line broadening in the chromosphere of a late type star. If an emerging loop were stable, then it would contribute little or nothing to macroturbulent velocity fields. Therefore, it is an essential part of the current picture that the magnetic flux loops in RS CVn stars and late type giants are unstable, moving upwards rapidly through the stellar chromosphere, with material draining down along both legs for at least some time during the evolution. The macroturbulent velocity parameter is, in this scenario, to be interpreted as related to the rise velocity of the unstable loops. In this regard, the sun provides a useful calibration: young active regions, where new flux loops are continually emerging, have been found to be the site of large macroturbulence (Shine and Linsky, 1973), and the value of the macroturbulence broadening parameter, ζ, in the brightest active regions ($\zeta \approx 10$ km/sec) agrees quantitatively with the upward velocity of rising flux loops (Bruzek and Durrant, 1977). Older active regions show much reduced ζ (Shine and Linsky, 1973), and in these, loops are no longer emerging.

In our scenario, the interpretation of line broadening in a stellar profile in terms of macroturbulent broadening requires a large number of macroscopically large elements on the surface of the star. Recently, various authors have made estimates of the numbers of loops which are present in the visible atmosphere of RS CVn stars, using a scaling relation which was originally derived for static loops in the solar corona (Walter et al., 1980; Swank et al., 1981). These authors have found that the numbers of loops

required on the visible hemisphere to explain the observed X-ray emission from RS CVn stars are large, 10^3-10^4. Now, application of static loop scaling relations cannot be quantitatively accurate if our scenario of non-static loops holds true. Thus, we cannot interpret these loop numbers as literally accurate. However, they are large enough to suggest that even if the condition of static loops were relaxed, the numbers of loops would still be large.

We are therefore led to propose that the macroturbulent "elements" which give rise to the large line broadening in λ And and other late-type stars are to be identified with unstable magnetic flux loops. Each loop would contain an essentially distinct "chromosphere", shielded by the magnetic field lines from the ambient medium through which it was passing. Hence, there is no difficulty with the fact that the macroturbulence velocities which we have been discussing (30-40 km/sec) are highly supersonic in the chromosphere: the magnetic field will shield the "chromosphere" in the rising loop from the effects of supersonic turbulent dissipation in the ambient medium.

If macroturbulence is determined by flux loops, it follows that the broadening parameter, ζ, is intrinsically time-dependent, with a value which depends on the immediate pre-history of flux loop emergence prior to the time of observation. At times, the number of loops will be small, in which case, it is expected that a line profile might break up into separate components, one from each loop. In fact, the Ca K line profile shows this kind of behavior (Baliunas and Dupree, 1979): at times it is flat topped (implying large numbers of rapidly moving macroturbulence elements), and at other times, a strong central absorption develops, indicating that the macroturbulence broadening has dropped to a small value.

The existence of a finite number of loops on the surface also results in some asymmetry in the line profiles. From purely statistical arguments, we can estimate that with 10^3-10^4 loops on the visible disk at any one time, the variations due to random changes in numbers of flux loops are expected to be of order 1-3%. Asymmetries of just this order are known to occur both in Ca K emission and in the Hα line profile in λ And (Baliunas and Dupree, 1979; Mullan and Cram, 1982).

(viii) The magnetic key: RS CVn stars

Our proposal for mass loss in cool giants rests on the assumption that there are magnetic fields in these stars. Nothing in the literature, however, forces us to believe unambiguously that late-type giants and supergiants have magnetic fields on the magnitude required to drive rapid mass loss in a series of discrete reconnection events. On the other hand, it seems rather likely that magnetic fields will exist in cool giants because of dynamo action in their convection zones. However, in order to put our proposal on a firm footing, we need to find a group of stars in which the proposed mechanism is at work, and which are known from other reasons to possess magnetic fields in their atmospheres. RS CVn active secondaries satisfy this requirement: circular polarization in their radio emission, and large starspots on their surfaces, provide strong circumstantial evidence for magnetic fields. The various pieces of evidence which have been accumulated in many spectral regions for these systems suggest flaring activity is in progress. Estimates of current rates of mass loss (Walter et al, 1978; DeCampli and Baliunas, 1979) seems at first sight to be small (few times 10^{-9} M_{sun}/yr). However, it is important to

realize that the average rate of mass loss and/or exchange between components over evolutionary time scales must be much smaller than this (10^{-11} M_{sun}/yr; Popper and Ulrich, 1977). Hence, the current mass loss process cannot have been in existence for more than about 1% of the stellar lifetime: i.e. it is a very recent development. This is consistent with our discovery that active secondary stars in RS CVn systems lie very close to the mass-loss boundary (as seen in Mg II h and k; Mullan, 1981).

Thus, all of the ingredients (flaring activity, recent mass loss, magnetic fields) appear to be present in the RS CVn stars, and we therefore consider these systems as the key to understanding the relationship between rapid mass loss in cool giants and magnetic activity. The RS CVn systems even provide evidence for discrete ejections of mass (Pfeiffer, 1979).

The reason that the RS CVn stars stand out against the background of all other late-type giant stars is that they have magnetic fields which are strong enough to make magnetic activity detectable at Earth. (Membership in a binary probably helps to make the magnetic fields strong by means of tidal action; Mullan, 1975). We propose that other late-type giant stars also have magnetic fields in their atmosphere, not strong enough in general to produce magnetic activity at a high enough level to be detectable readily at earth, but still strong enough to have a dominant dynamic effect on mass loss from that atmosphere. And although the flaring activity in these other late-type stars is not as prominent as in RS CVn systems, nevertheless, evidence is beginning to accumulate which shows that some activity is indeed in progress in late-type giants (cf. Wischnewski and Wendker, 1981; Boice et al., 1981; Baliunas et al., 1981; Mullan and Stencel, 1982).

(ix) C^{13} richness

Perhaps the most interesting feature to emerge in the discussion of "boundary lines" in the HR diagram in recent months involves the heavy isotope of carbon, C^{13}. This isotope is created by CNO burning inside massive stars. During evolution of these objects, it must presumably be dredged up from deep within the star if it is ever to appear in the atmosphere in amounts which exceed the "normal" abundance (i.e. the interstellar medium abundance of about 1/89 of C^{12}). There are severe problems with most of the mechanisms which have been devised so far to explain the required mixing, not only of C isotopes, but also of other elements (cf. Iben, 1981; Scalo, 1981). After a long discussion of many possible mixing scenarios, Scalo (1981) concludes "our understanding of the advanced stages of stellar evolution is seriously incomplete, a stochastic mixing process is at work" (Scalo, 1981, p. 104).

It is with this in mind that we draw attention to an important feature of C^{13} rich stars: <u>they show a strong preponderance to lie on the rapid mass loss side of the Mg boundary</u> discovered by Stencel and Mullan (1980). Thus, in a sample of stars studied by Lambert and Ries (1981) for C^{13} abundance, 18 stars overlap with our sample of mass loss stars. Of these, 11 show S/L > 1 in our data (no rapid mass loss), and 7 show S/L < 1. Of the former, 9 have S/L > 1 in our Mg data (i.e. solar-like atmospheres, no rapid mass loss), while of the latter, 5 have S/L < 1, while the other 2 have S/L $\approx$ 1. Thus, <u>C^{13} richness appears to be rather highly correlated with rapid mass loss</u>.

How can this be fitted into our scenario of unstable magnetic loops? Obviously, the unstable loops are dredging up material from deep within the star, wherever the dynamo action is occurring. It is believed (Rosner, 1980) that dynamo activity occurs rather deep inside many stars, namely, beneath the convection zone. In red giants, this would place the dynamo region certainly deep enough to have access to CNO-processed material. In fact, magnetic mixing has already been proposed as an explanation for elemental composition in one of the stars where mass loss is rapid (Arcturus: Hubbard and Dearborn, 1980), and Scalo (1981) mentions magnetic fields as a possible candidate for mixing of elements in peculiar giants in general. The characteristic of stochasticity required for the mixing processes (Scalo; quoted above) is obviously well satisfied by a magnetic loop mechanism, since loops are known to appear at the surface in a chaotic fashion in time.

IV. T TAURI STARS

I have used empirical estimates of pressures in T Tau chromospheres (Cram, private communication) to estimate that these stars lie close to the MTTL, just as active secondaries of RS CVn stars do. However, there is one important distinction: RS CVn stars are evolving across the MTTL from left to right (they are older stars, 10^9 years old, on the basis of tidal synchronization; Mullan, 1981) while T Tau stars are young and are evolving in the opposite direction. Thus, there seems to be no difficulty in ascribing high macroturbulence, activity and mass outflow rates in T Tau stars to the same physical mechanism as in RS CVn stars i.e. unstable magnetic loops.

The distinction about the direction of evolution, however,

makes for some differences between RS CVn stars and T Tau stars. For example, C^{13} richness ought not to be a characteristic of rapid mass loss in T Tau stars, according to our scenario, since, evolution has not yet been able to proceed through CNO burning in T Tau stars. Moreover, since the T Tau stars are moving from right to left across the MTTL, they are moving in the direction of having magnetic loops become stable (rather than unstable). Hence, we expect that some loops, rather than breaking away from the star, may evolve outwards for a time, and then be restrained by the stellar gravity, and become a stable loop. Then, upward motion would cease, and only material draining back down along the loop legs would be visible in the spectrum, and material would be observed to be falling into the star. Thus, T Tau stars ought to be characterized by both outflow and inflow of material, depending on whether the loop is on the unstable side of MTTL, or on the stable side. If most T Tau stars are on average above MTTL, then the lack of correlation between chromospheric and coronal emission reported by several authors (e.g. Walter and Kuhi, 1981) can be understood in terms of the arguments given above (section III (iii)).

V. CONCLUSION

Unstable emerging magnetic flux loops have emerged as a unifying factor in our study of RS CVn systems. However, although the idea is most strongly supported among the RS CVn systems, the idea appears to have much broader application, namely, it may be applicable to most late-type giants which are losing mass rapidly. We have been able to use the concept of unstable loops to account for rapid mass loss, episodic mass loss, flaring activity, high macroturbulence in chromospheres, He 10830 A line data (i.e. maximum

strength correlated with maximum activity), and C^{13} richness in rapid mass loss stars.

The work which has been done in the course of this grant has opened up several avenues for future research, avenues which have an impact on the way in which we will approach not only the question of mass-loss and activity, but also the "solar-stellar connection" in general.

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FIGURE CAPTIONS

Figure 1. Examples of Hα flux profiles from various stellar model chromospheres. The photospheric model is specified in the lower left corner. G denotes that it is model of Gingerich (Carbon and Gingerich, 1969), the first three numbers denote $T_{eff}/10$, and the last three are 100 log g. The four curves for each photospheric model refer to four chromospheres, with T_c, log m_c values as follows: _._._._ 6500, -6; _.._.._.._.. 6500, -4; ----- 7500, -4; .... 7500, -6. In the model with T_{eff} =6000 K, log g = 2.00, the peculiar behavior in one curve is due to data drop out during plotting.

Figure 2. Observed and theoretical line profiles of Hα in λ And. Observations refer to 2 of 4 epochs provided to us by Bopp and Smith (1980). Model 1 is our low pressure chromosphere, broadening by macroturbulence; Model 2 is a high pressure chromosphere, without turbulence. The observed profiles have been extracted from the data of Bopp and Smith within 2Å of line center: the observed line is somewhat asymmetric, and is not completely smooth; we have chosen the smoother side of line center to extract the half profiles shown here.

Figure 3. Scenario for magnetically driven mass loss. A new flux loop emerges at time t_0 from beneath the stellar surface, and balloons upwards. Reconnection near the base eventually leads to a severing of the magnetic link with

the star, and a bubble of material is ejected. Each new flux loop which becomes unstable contributes a new mass loss episode in this scenario.

Figure 4. Survey of chromospheric velocity fields in the HR diagram shows a dividing line (VDL=velocity dividing line) between stars where mass loss is rapid (X) from those where mass loss rates are small (o) (Data from Stencel and Mullan, 1980). TDL denotes temperature dividing line (Linsky and Haisch, 1979; see also Simon et al., 1982 for further confirmation of the existence of TDL.) Note that cool coronae set in at essentially the same place where mass loss becomes rapid. Circled P's and π's denote STL (supersonic transition locus; Mullan, 1978) and MTTL (magnetic topology transition locus; Mullan, 1981, 1982) according to evolutionary tracks of Paczynski (1970). Note that MTTL and VDL are essentially coincident, i.e. mass loss becomes rapid in those stars where static closed loops of magnetic flux can no longer exist.

PERSONNEL

During the period of this grant, Dr. Wayne Waldron was appointed as a post-doctoral fellow at Bartol. Dr. Waldron has a master's degree in Plasma physics from Ohio State University, a master's degree in plasma physics from University of Wisconsin, and a Ph.D. in Astrophysics from the University of Wisconsin. His Ph.D. work was on winds from hot stars, the effects of hot coronae, and how these affect the radiatively driven winds in these stars. His work has involved an important new step in analyzing X-rays from hot stars, for his model allows one to follow the transfer of X-rays outwards from a base corona through a wind. Thus, he has been able to include an energy equation to follow how the effects of recombination and photo-ionization control the temperature structure of the wind. His detailed knowledge of stellar winds, and also of MHD effects in plasma physics, make him an ideal person to undertake the problem described in this report on magnetically driven mass loss from stellar coronae. He has been working to develop a 2-D MHD code to study the ballooning instability of a flux loop emerging into a stratified atmosphere.

PAPERS PRESENTED AT MEETINGS:

1. "Non-thermal Stellar Winds in Cool Stars", presented by D. J. Mullan at Cool Star Workshop, Cambridge, MA, January 31, 1980.
2. "Mass Loss from Warm Giants: Magnetic Effects", presented by D. J. Mullan, Erice Workshop on Physical Processes in Red Giants, September 3-13, 1980.
3. "Variable Mass Loss and Magnetic Topology in Cool Giant Stars", by D. J. Mullan and R. E. Stencel, presented by D. J. Mullan, American Astronomical Society Meeting, Calgary, June-July 1981.

4. "Large Macroturbulence in the Chromosphere of an RS CVn Star", by L. E. Cram and D. J. Mullan, presented by D. J. Mullan, American Astronomical Society Meeting, Calgary, June-July, 1981.
5. "Are discrepant asymmetry red giants necessarily hybrid stars?" D. J. Mullan and R. E. Stencel, presented at American Astronomical Society Meeting, Boulder, January 1982.
6. "Discrepant Asymmetry Stars: the Role of Unsteady Magnetic Flux Loops in the Atmospheres of Late-Type Giant Stars", D. J. Mullan and R. E. Stencel, to be presented at "Four Years of IUE", March, 1982.
7. "Static and Non-Static Helmet Streamers in Stellar Atmospheres: Effects on Mass Loss in Cool Stars", D. J. Mullan, to be presented at IAU Symposium No. 102, "Solar and Stellar Magnetic Fields: Origins and Coronal Effects", Zurich, August 1982.

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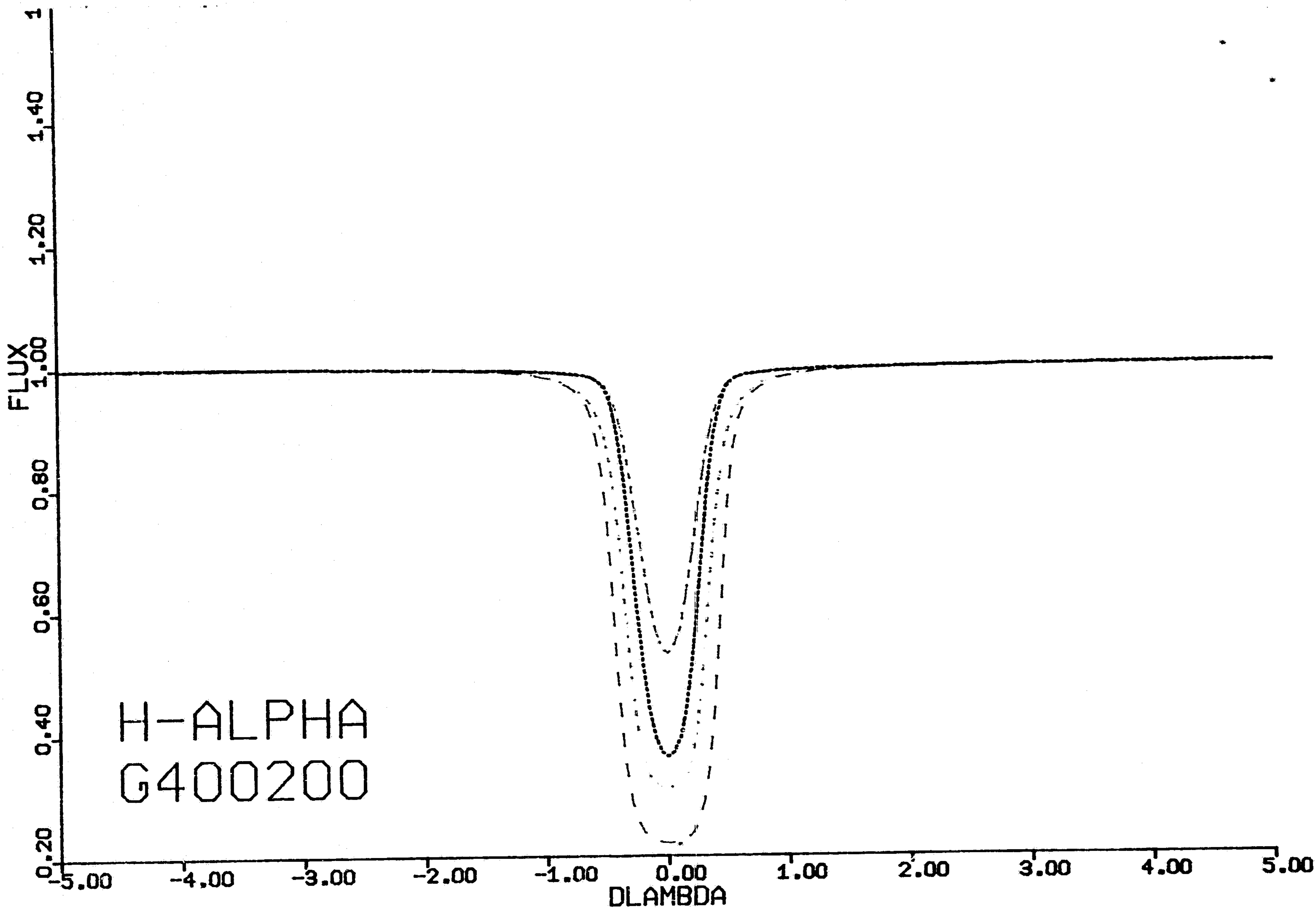
FLUX
1
1.40
1.20
1.00
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-5.00
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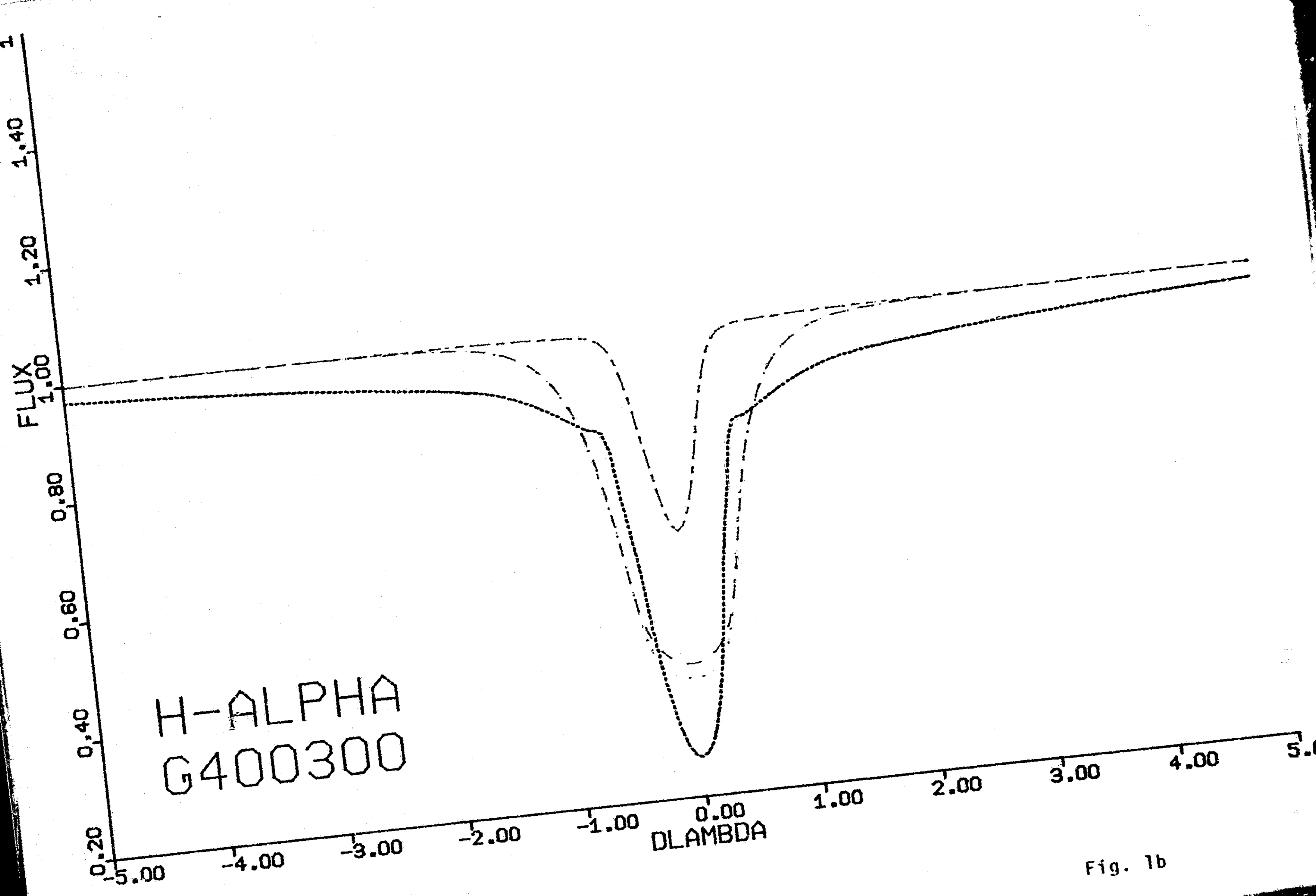
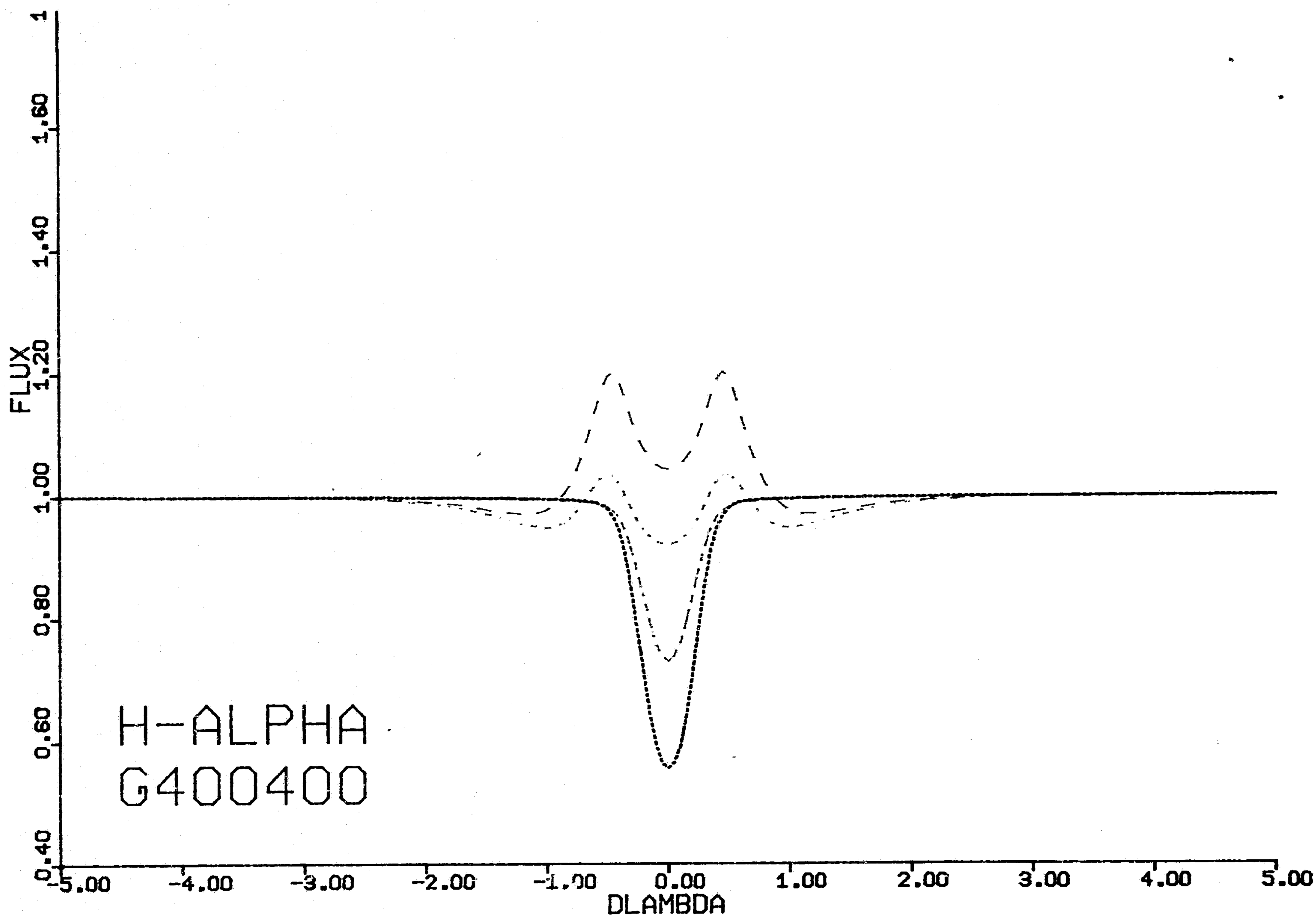
G400200

Fig. 1a



H-ALPHA
G400300
FLUX
1
1.40
1.20
1.00
0.80
0.60
0.40
0.20
DLAMBDA
-5.00
-4.00
-3.00
-2.00
-1.00
0.00
1.00
2.00
3.00
4.00
5.00

Fig. 1b


H-ALPHA
G400400
FLUX
DLAMBDA
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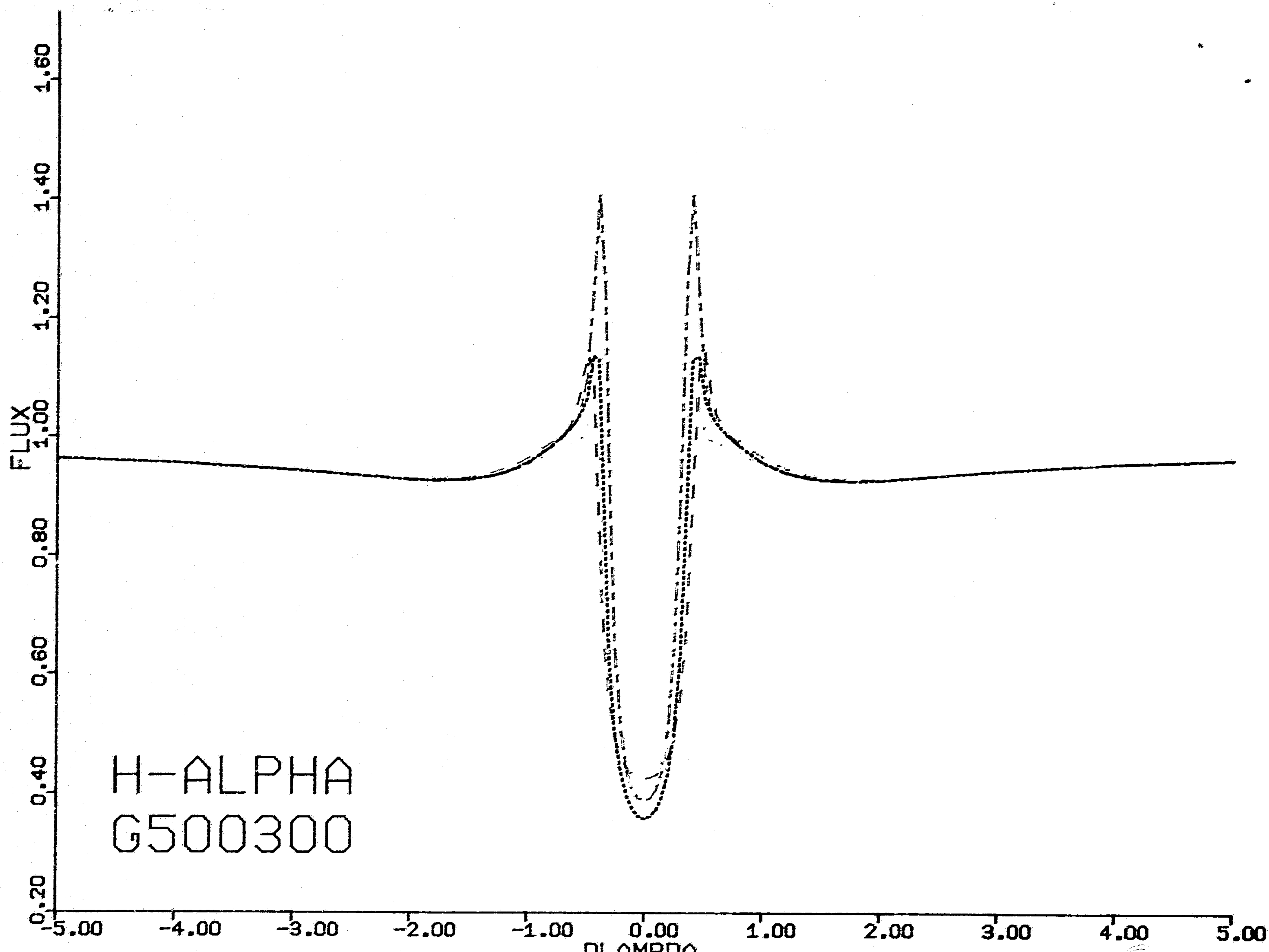
0.60
0.80
1.00
1.20
1.40
1.60
1
-5.00
-4.00
-3.00
-2.00
-1.00
0.00
1.00
2.00
3.00
4.00
5.00

Fig. 1c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G500300
FLUX
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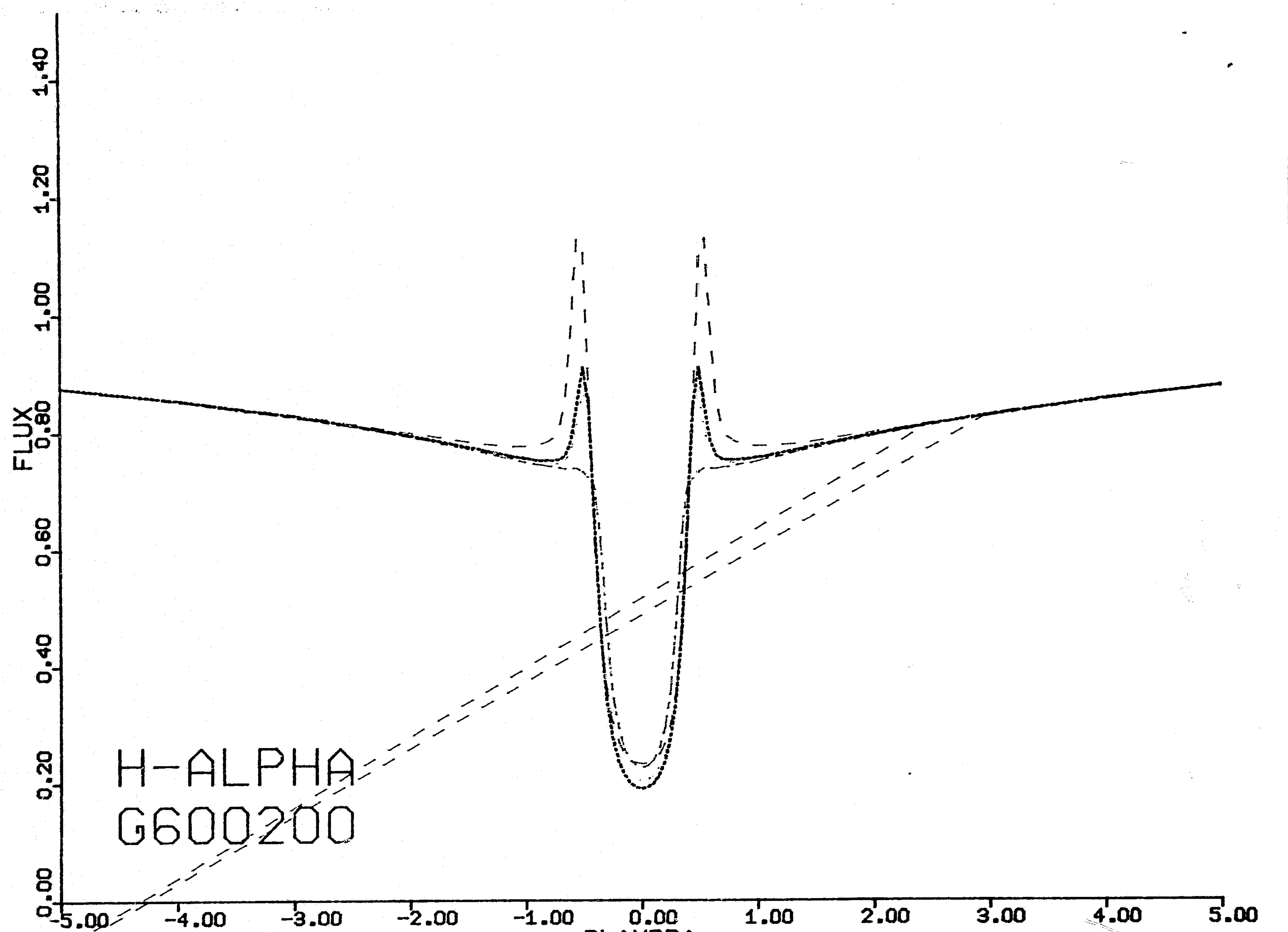
1.40
1.20
1.00
0.80
0.60
0.40
0.20
-5.00
-4.00
-3.00
-2.00
-1.00
0.00
1.00
2.00
3.00
4.00
5.00


H-ALPHA
G600200
FLUX
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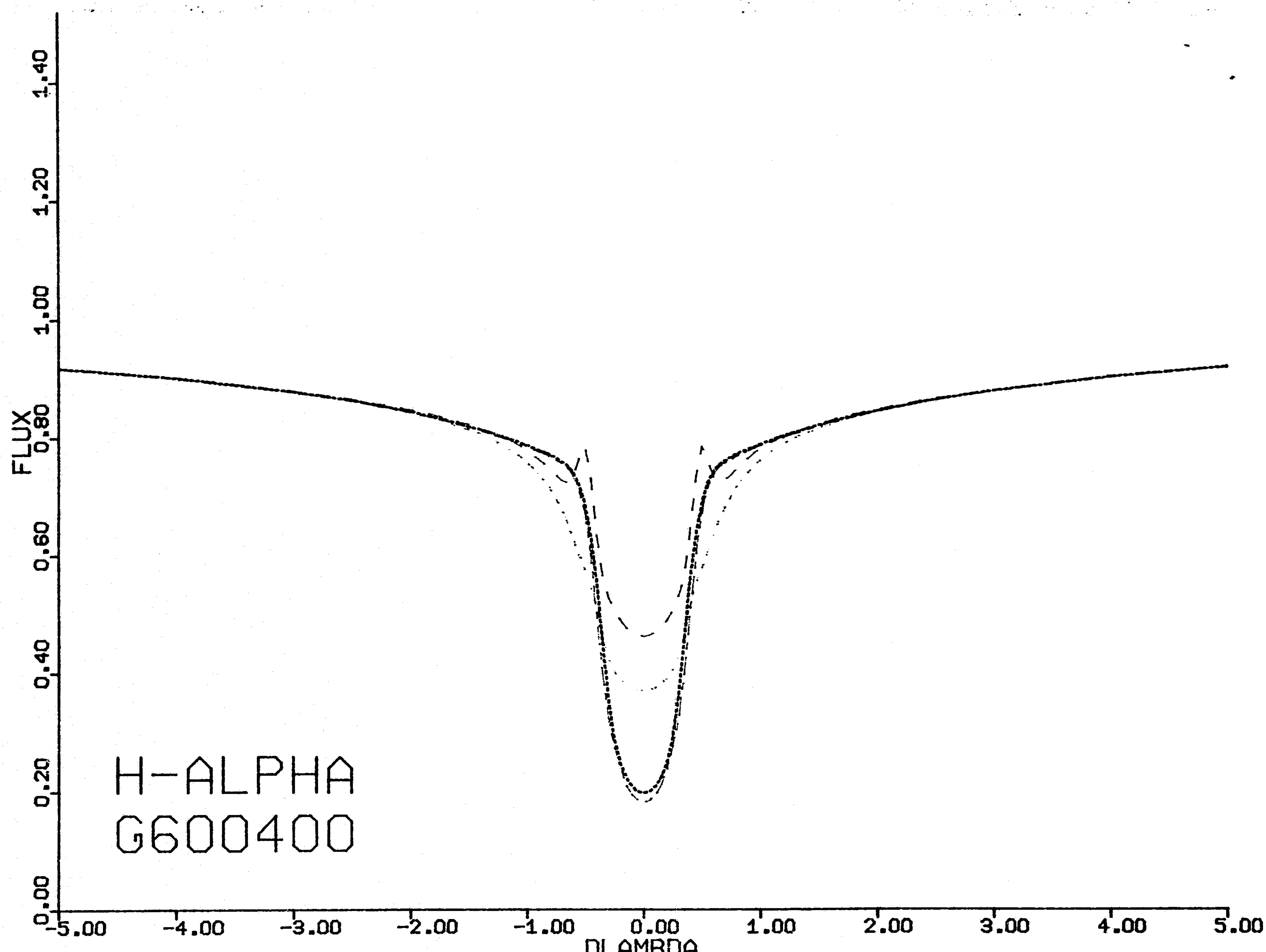
1.20
1.00
0.80
0.60
0.40
0.20
0.00
-5.00
-4.00
-3.00
-2.00
-1.00
0.00
1.00
2.00
3.00
4.00
5.00



H-ALPHA
G600400
FLUX
1.40
1.20
1.00
0.80
0.60
0.40
0.20
0.00
-5.00
-4.00
-3.00
-2.00
-1.00
0.00
1.00
2.00
3.00
4.00
5.00
DLAMBDA

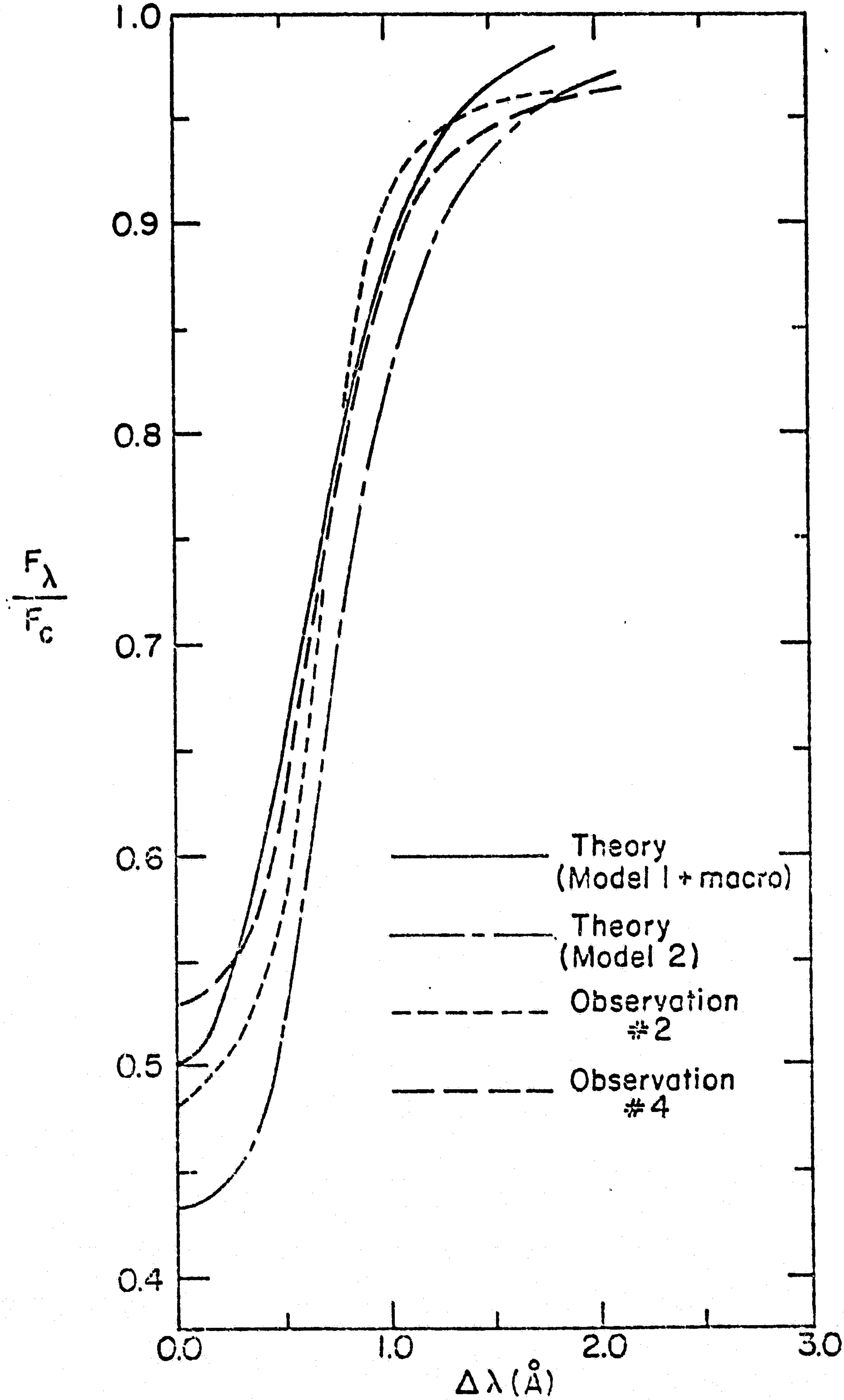

1.0
0.9
0.8
0.7
0.6
0.5
0.4
$\frac{F_\lambda}{F_c}$
0.0
1.0
2.0
3.0
$\Delta\lambda$ (Å)
Theory (Model 1 + macro)
Theory (Model 2)
Observation #2
Observation #4

Fig. 2

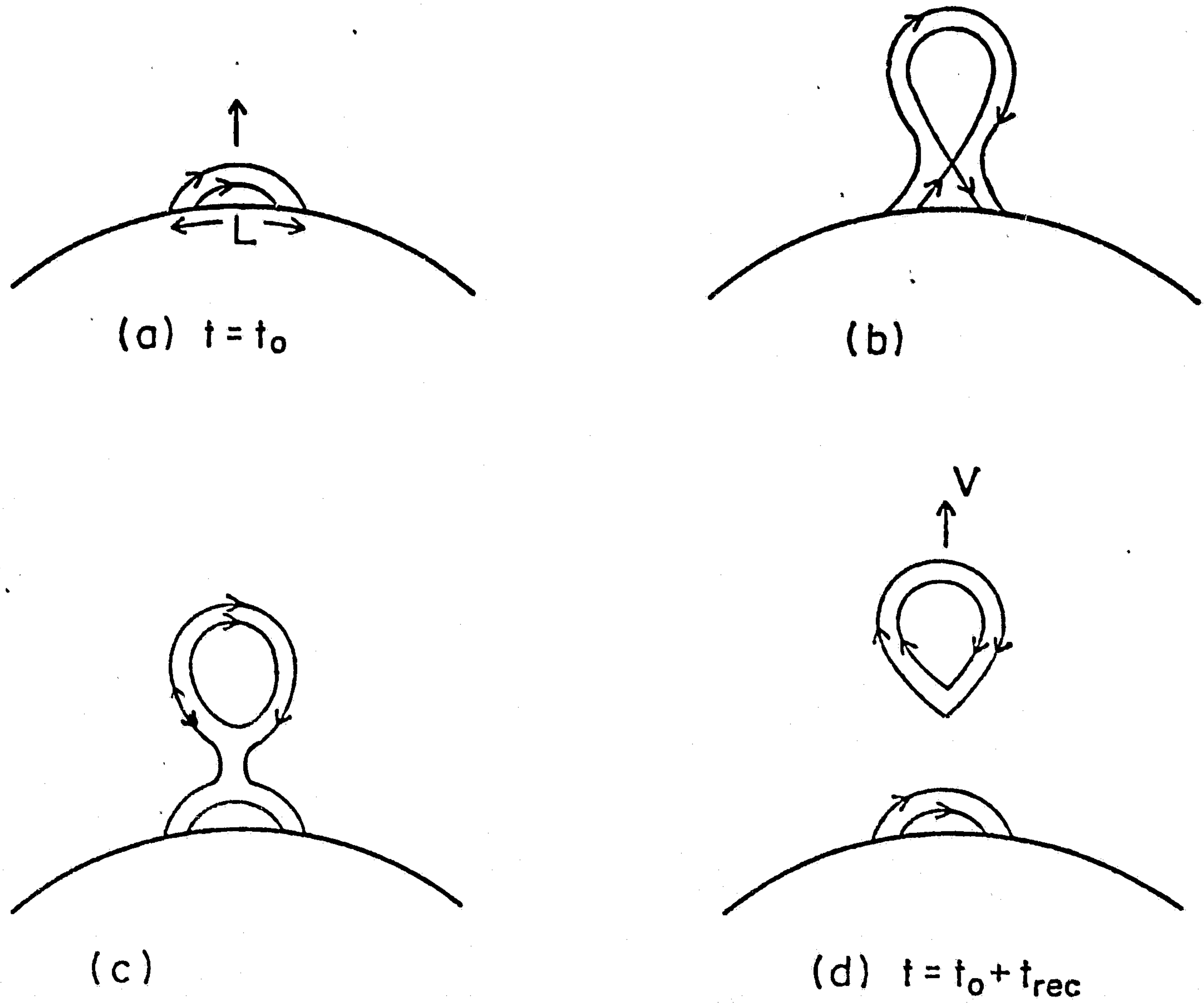


Fig. 3

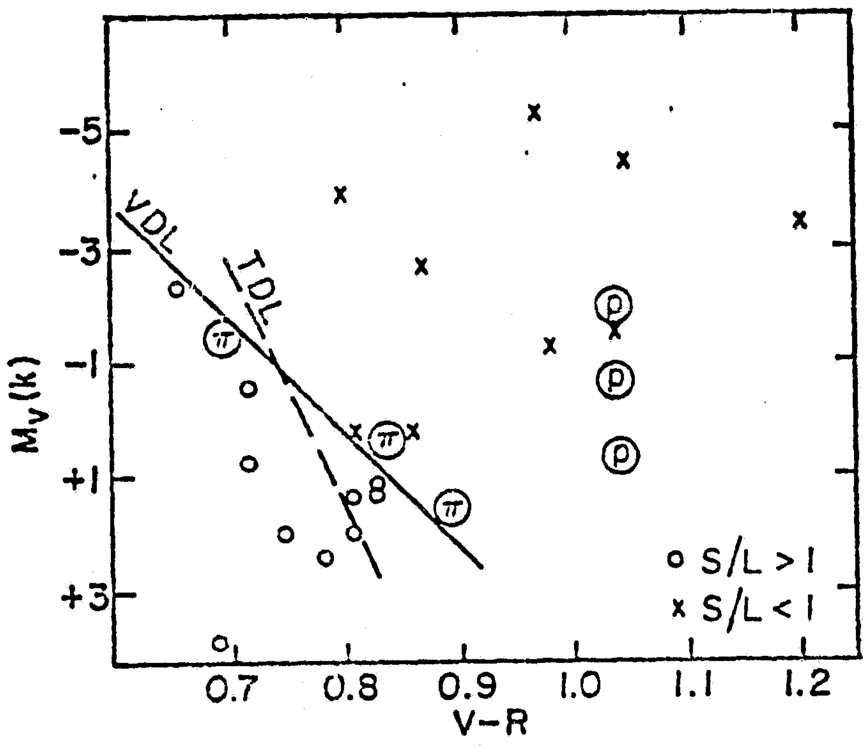


Fig. 4